B[E] SUPERGIANTS: WHAT IS THEIR EVOLUTIONARY STATUS?

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Abstract. In this paper, we investigate the evolutionary status of B[e] stars from the point of view of stellar evolution theory. We try to answer to the question of how massive hot supergiants — i.e. evolved stars — can be capable of producing a circumstellar disk. We find and discuss three possibilities: very massive evolved main sequence stars close to critical rotation due to their proximity to their Eddington-limit, blue supergiants which have just left the red supergiant branch, and single star merger remnants of a close binary system. While the latter process seems to be required to understand the properties of the spectroscopic binary R4 in the LMC, the other two scenarios may be capable of explaining the distribution of the B[e] stars in the HR diagram. The three scenarios make different predictions about the duration of the B[e] phase, the time integrated disk mass and the stellar properties during the B[e] phase, which may ultimately allow to distinguish them observationally.

1. Introduction

In recent years, the so called B[e] supergiants (cf. Lamers, this volume) have emerged as a distinct class of massive stars, whose defining properties — a strong mid-IR excess, strong Balmer emission, and narrow permitted and forbidden low-excitation emission lines — can be explained by a slowly outflowing equatorial disk superimposed to a normal fast wind (cf. Zickgraf et al. 1996a). Although the number of these objects is small — the best statistics exists for the Magellanic Clouds, where about 15 B[e] supergiants

have been found (Gummersbach et al. 1995) — it is comparable to the number of Luminous Blue Variables (LBVs; cf. Bohannan 1997), which are know to represent a key evolutionary phase of very massive stars (Langer et al. 1994, García-Segura et al. 1997, Langer et al. 1998). It is thus of fundamental importance to understand whether the B[e] supergiants are just freaks, i.e. peculiar objects which come to exist due to special circumstances — in which case they might still show interesting physical phenomena — or whether all stars within a certain initial mass range evolve through a B[e] phase. In this case, our general understanding of the evolution of stars in that mass range may depend on our understanding of the B[e] stars.

Although we have a solid knowledge of the evolution of the deep interior of massive stars since a long time (Weaver et al. 1978, Kippenhahn & Weigert, 1990) it has become more and more clear during the last decades that our understanding of the evolution of their observable features is still rather incomplete. Among others, open problems concern the effective temperature evolution of moderately massive ($M_{\rm ZAMS} \simeq 10...30\,M_\odot$) post main sequence stars (Langer & Maeder 1995) and the evolutionary connections between O stars, Luminous Blue Variables (LBVs) and Wolf-Rayet (WR) stars for higher masses (Schaller et al. 1992, Langer et al. 1994, Stothers & Chin 1996, Pasquali et al. 1997; see also Schulte-Ladbeck 1998). Two major physical difficulties in the theoretical models of the observable evolutionary stages of massive stars have been identified and made responsible for the persisting lack of reliable models: mass loss and internal mixing processes (Meynet et al. 1994, Langer 1994, Deng et al. 1996). E.g., it has been found that the mass loss of massive main sequence stars should be roughly twice as high as what appears to be observed in order to understand many features of massive post main sequence stars (Meynet et al. 1994, Langer et al. 1994). Additionally, there is growing evidence that stellar rotation may considerably affect the evolution of massive stars (Maeder 1987, Langer 1991a, Fliegner et al. 1996, Maeder & Meynet 1996, Meynet & Maeder 1997, Langer et al. 1997ab). *Rapid rotation* can reduce the effective gravity in the star, and it produces large scale flows (Eddington 1925). During the evolution, *differential rotation* occurs in all stars, with the possibility of the occurrence of various local hydrodynamic instabilities (cf. Endal & Sofia 1978, Zahn 1983) and corresponding mixing of chemical elements and angular momentum. Of relevance for massive stars are the shear instability (cf. Maeder 1997), the baroclinic instability (Zahn 1983, Spruit & Knobloch 1984), and the Solberg-Høiland and Goldreich-Schubert-Fricke instabilities (cf. Korycansky 1991).

Time dependent evolutionary models for massive stars including rotation have been constructed in the past in one dimension, using various degrees of approximation (e.g. Endal & Sofia 1978, Maeder 1987, Langer

1991a, Langer 1992, Talon et al. 1997, Langer 1998; cf. also Dupree 1995). Today, it is beyond reasonable doubts that the evolution of massive stars is influenced by rotation due to the physical mechanisms mentioned above (cf. Fliegner et al. 1996). While the principle effects of rotation in the interior of massive stars during their evolution all the way to iron core collapse are described elsewhere (Langer et al. 1997b, Heger et al. 1998), we discuss here whether massive single stars can approach the limit of critical rotation. We shall assume in Sections 2 and 3 that this would lead to the disk structure responsible for the B[e] phenomenon, either due to the Bjorkman-Cassinelli (1993) mechanism of wind compression of a rotating star or otherwise. A detailed wind disk model for B[e] stars can be found in Bjorkman (this volume), while the recent status of the Bjorkman-Cassinelli model is discussed by Owocki et al. (1996), Owocki (this volume) and Cassinelli (this volume). In Section 4 we will discuss a model where the disk is not produced by a critically rotating single star but rather a critically rotating close binary.

2. Very massive main sequence stars

Massive main sequence stars are rapid rotators, with equatorial rotation velocities in the range of 100...400 $\mathrm{km\,s^{-1}}$ (Fukuda 1982, Penny 1996, Howarth et al. 1997). In our first scenario, we investigate whether massive main sequence stars, i.e. massive stars during core hydrogen burning, are capable of arriving a critical rotation (cf. Langer 1998). The results of this Section have been obtained with a hydrodynamic stellar evolution code (cf. Langer et al. 1988, Langer 1991b), employing the OPAL opacities of Iglesias et al. (1992). We have used an outer boundary condition which takes the optical depth of the stellar wind selfconsistently into account within a grey approximation (Langer et al. 1994, Heger & Langer 1996).

We have computed the proximity of the star to the Eddington-limit, i.e. the Eddington factor $\Gamma = L/L_{\rm edd} = \kappa L/(4\pi c G M)$ in the following way. It has been shown in Langer (1997) that the occurrence of convection and of density inversions makes the concept of the Eddington limit as a stability limit invalid in the stellar interior, i.e. the Eddington factor Γ has to be evaluated only at the stellar surface. Since the term “surface” is not unambiguously defined in this context, we considered Γ in layers with an optical depth of $\tau < 100$, where in fact neither a significant convective energy flux nor density inversions have been found in the investigated models. Thus, to estimate the distance to the Eddington limit, we used the maximum value of Γ occurring in the subsonic layers with $\tau < 100$. Furthermore, we used the OPAL opacity coefficient to compute the Eddington factor. Mass and luminosity are practically constant for $\tau < 100$ and equal to the total stellar mass and luminosity.

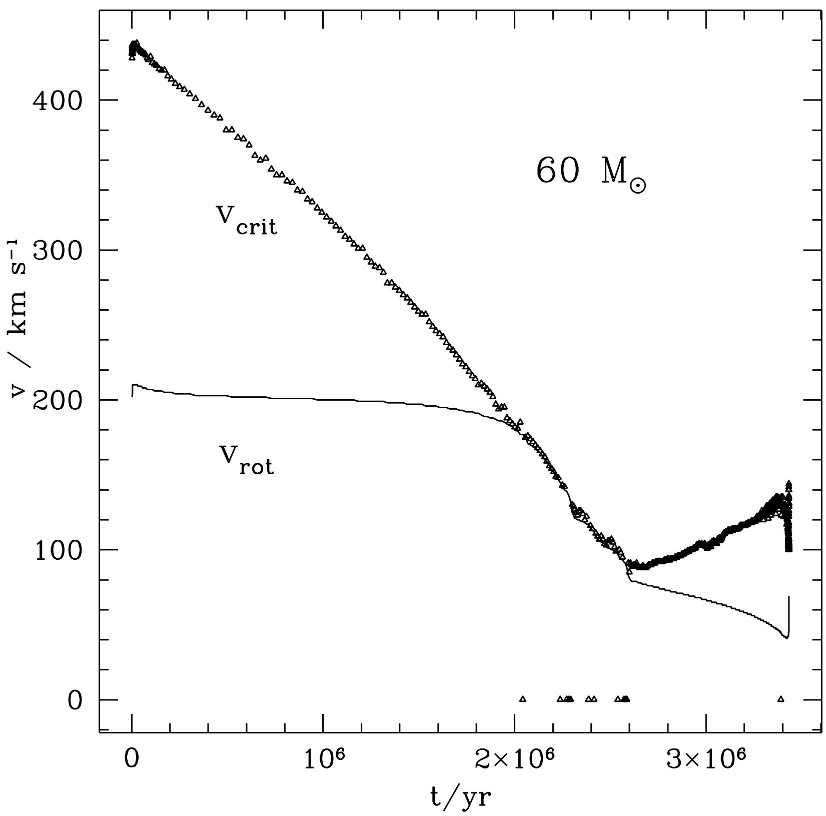

Figure 1. Equatorial rotational velocity $v_{\rm rot}$ and the critical rotational velocity $v_{\rm crit}$ as function of time during the main sequence evolution of a 60 $M_\odot$ sequence.

Angular momentum is carried only as a passive quantity in the stellar models, i.e. we ignore the centrifugal force in the stellar interior as well as the effect of mixing of chemical species due to rotationally induced instabilities (cf. Fliegner et al. 1996, Meynet & Maeder 1997). However, we do consider the centrifugal force at the stellar surface to evaluate the distance of the star from the Ω-limit (Langer 1997), i.e. from critical rotation, with $\Omega = v_{\rm rot}/v_{\rm crit}$, and $v_{\rm crit}^2 = GM(1-\Gamma)/R$. Furthermore, we assume our models to be always rigidly rotating. Stellar models including differential rotation (e.g. Fliegner et al. 1996) show that this is a good approximation on the main sequence, since the time scale for angular momentum transport is of the order of the thermal time scale and also shorter than the

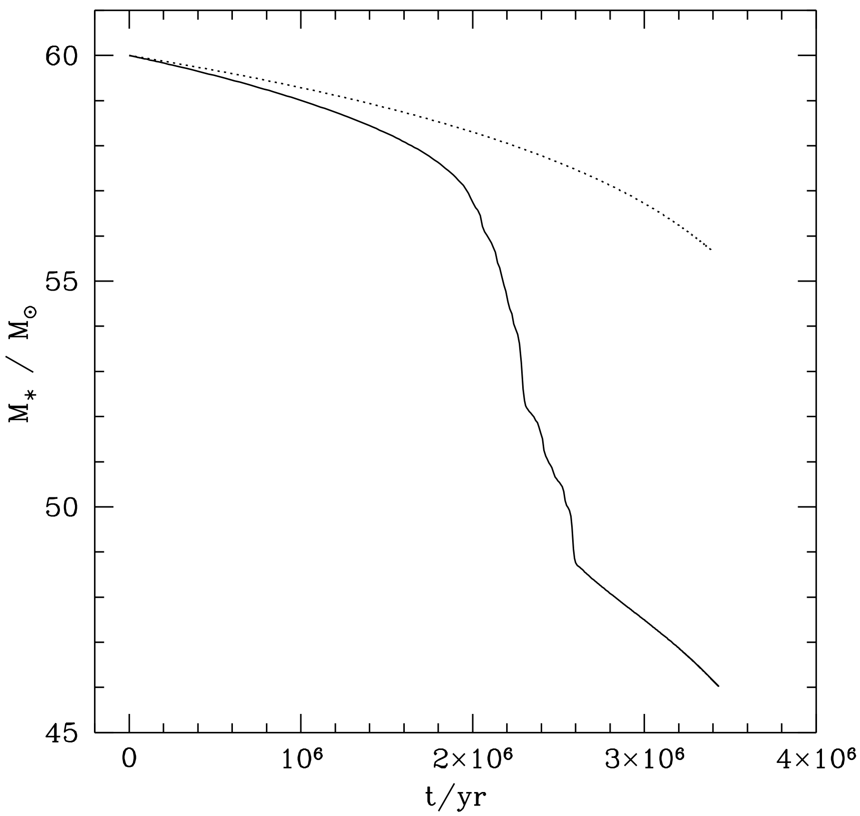

Figure 2. Stellar mass as a function of time for the $60\,M_{\odot}$ sequence shown in Fig. 1 (solid line), in comparison to the evolution of a $60\,M_{\odot}$ star without enhancement of the mass loss rate due to rotation (dotted line). The mass loss rates can be read of this figure as the slope of the curves.

time scale of rotationally induced chemical mixing (Chaboyer & Zahn 1992, Zahn 1992, Talon & Zahn 1997). According to Zahn (1994), the expected amount of differential rotation in a massive main sequence star is roughly $\Delta\omega/\omega \simeq \omega^2 R^3/(GM)$, with ω being the mean angular velocity and $\Delta\omega$ its difference between stellar core and surface. In the models presented below, this estimate gives $\Delta\omega/\omega \lesssim 0.01$. The approximation of rigid rotation may become invalid when the mass loss time scale becomes shorter than that of angular momentum transport, but this is not the case in our models.

Angular momentum loss is only considered through the effect of mass loss, i.e. the lost mass carries away its specific angular momentum. To com-

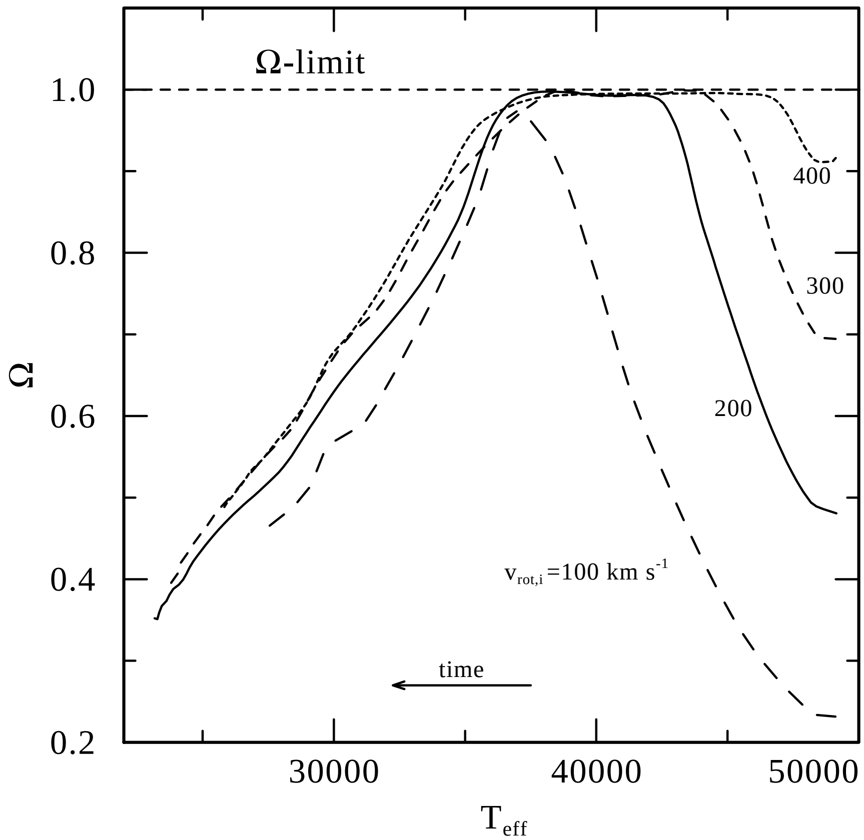

Figure 3. Evolution of the ratio of the rotation rate to the critical rotation rate, Ω, as function of the effective temperature during core hydrogen burning, for 60 $M_\odot$ sequences starting with different equatorial rotational velocities $v_{\rm rot,i}$.

pute the mass loss rate for our stellar models, we have applied the empirical rate found by Lamers & Leitherer (1993) with the metallicity dependence obtained by Leitherer & Langer (1991). However, we have applied the correction factor derived by Bjorkman & Cassinelli (1993) as a fit to the results of Friend & Abbott (1986) to take the effect of rotation on the mass loss rate of hot star winds into account (cf. Langer 1998).

Fig. 1 shows the time evolution of the critical rotational velocity and the actual rotational velocity for a 60 $M_\odot$ sequence starting with $v_{\rm rot} = 200\,{\rm km\,s^{-1}}$. The critical rotational velocity has a pronounced minimum at roughly $t = 2.6\,10^6$ yr, which corresponds to a stellar effective temperature of $T_{\rm eff} \simeq 36\,500\,$K, around which the iron opacity peak has its maximum effect. We see in Fig. 1 that, as $v_{\rm rot}$ approaches $v_{\rm crit}$ (i.e. $\Omega \to 1$) the rotation

rate of the star declines such that the Ω-limit $\Omega = 1$ is never exceeded.

The reason is that, using the results of Friend & Abbott (1986), the mass loss rate increases as the star approaches $\Omega = 1$. In Section 3, we discuss that the angular momentum loss rate is directly coupled to the mass loss rate of the star (cf. Langer 1998, and Fig. 6 below). Consequently, the mass loss rate at the Ω-limit is determined by the angular momentum loss rate which is required to prevent the star from reaching $\Omega = 1$.

The result is that the star evolves along the Ω-limit until the critical rotation rate increases again, due to changes in the photospheric parameters (i.e., the opacity). For the considered example, the time spent at the Ω-limit is roughly $6\,10^5$ yr. Fig. 2 shows that during this time the mass loss rate is increased to roughly $10^{-5}\ M_{\odot}\ \mathrm{yr}^{-1}$, which is a factor of ~ 10 above the normal radiation driven mass loss rate (cf. Fig. 2).

According to this picture, very massive main sequence stars may thus, due to their proximity to their Eddington-limit, evolve at critical rotation for several 10^5 yr. Fig. 3 shows that, for a given star, this time scale depends on its initial rotation rate. As during the time at the Ω-limit the star may have a slow dense equatorial outflow, a relationship to the B[e] stars may be suggested.

3. Supergiants on a blue loop

In this Section we describe a possibility for a star to arrive at critical rotation which does not require an extremely high stellar luminosity. It was found to occur in contracting stars, whose envelope structure changes from convective to radiative (see Heger & Langer 1998). This situation occurs for massive stars undergoing a so called blue loop, i.e. core helium burning red supergiants which evolve off the Hayashi-line towards the regime of blue supergiants in the HR diagram. Blue loops are typically found in the initial mass range 5...25 $M_{\odot}$ (cf. Langer 1991b, Schaller et al. 1992). However, the mechanism discussed in the following may also apply to other situations, e.g. to the post-AGB phase of low mass stars.

Fig. 4 shows evolutionary tracks for rotating stars between 10 and 20 $M_{\odot}$ (cf. Heger et al. 1997 for details), two of which evolve through a blue loop. The 12 $M_{\odot}$ sequence will be investigated in more detail below. Figure 5 shows the time evolution of the radii of various Lagrangian mass shells and of the stellar radius, before and during the blue loop. Also shown is the radial extent of the convective part of the envelope. We see from Fig. 5 that, within a fair approximation, the inner and the outer radius of the convective envelope remain constant until the blue loop occurs, while mass shells continuously flow out of the convection zone through its lower boundary.

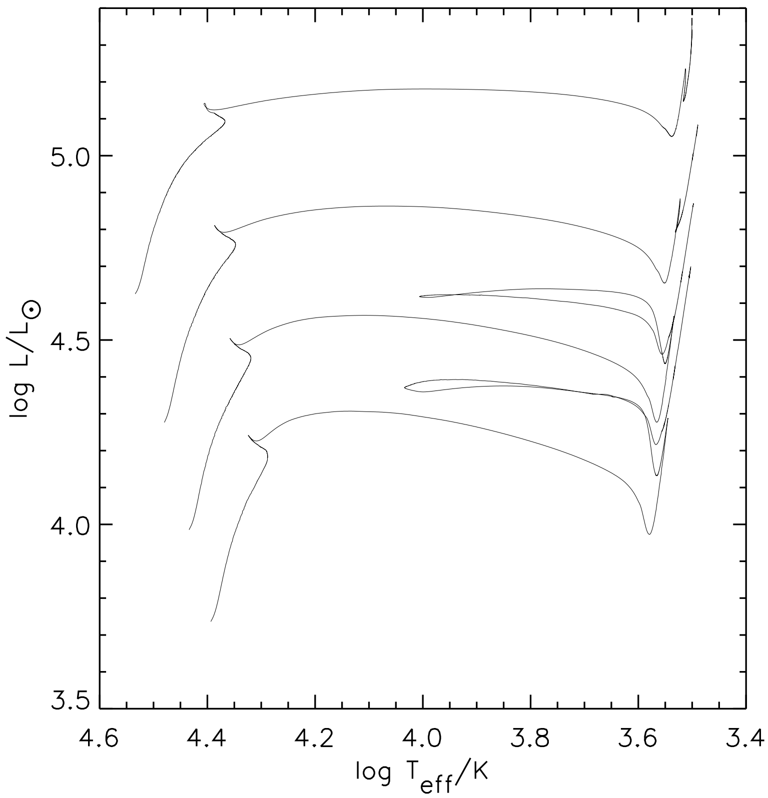

Figure 4. Evolutionary tracks for rotating 10, 12, 15 and 20 $M_\odot$ sequences from the ZAMS to central neon exhaustion (cf. Heger et al. 1997).

In Fig. 6 we show that, due to the rapid angular momentum transport within convection zones, this situation leads to a dramatic spin-up of the envelope. In the right part of Fig. 6 (B1...B4), the situation is sketched for the assumption of rigid rotation in convective regions and broken up into three discrete steps. For a constant inner radius of the convection zone, the drop-out of mass through its lower boundary leads to an increase of the rotation frequency and of the mean specific angular momentum in the convection zone. Note that rigid rotation is assumed only for simplicity; it is not necessary for the spin-up mechanism to work.

The left side of Fig. 6 (A1...A4) sketches the situation of the loss of mass through the upper boundary of a rigidly rotating region. In this case, which applies to the mass loss of very massive main sequence stars (cf. Sect. 2), matter with the highest specific angular momentum is continously

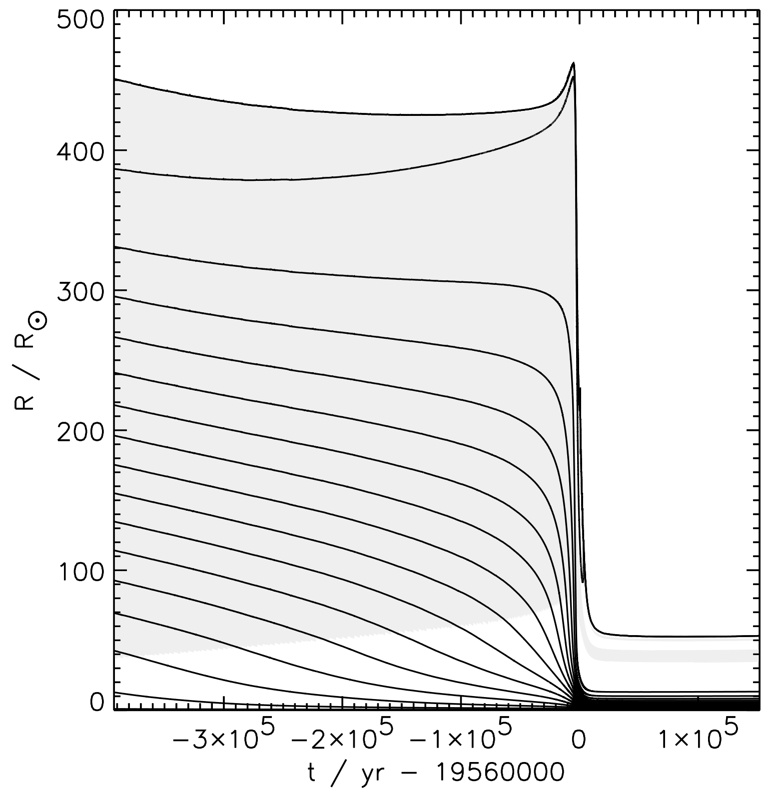

Figure 5. Evolution of the radii of different mass shells as a function of time for a period including the transition from the red to the blue supergiant stage of our 12 $M_\odot$ model. $t = 0$ is defined as in Fig. 7 and corresponds roughly to the time of the red-blue transition. Except for the uppermost solid line, which corresponds to the surface of the star, the lines trace Lagrangian mass coordinates. The mass difference between the lines is 0.5 $M_\odot$. Shading indicates convective regions.

lost, which leads to a spin-down of the rigidly rotating star.

Fig. 7 shows the time dependence of the rotational velocity of the 12 $M_\odot$ sequence before, during, and after the red-blue transition. The transition itself takes only about 10 000 yr. During this time the angular momentum transport to the surface layers of the star continues. The envelope layers remain rigidly rotating, continuing to shovel up angular momentum to the surface. The contraction of the star by a factor of $f \approx 10$ would increase the rotational velocity by the same factor if j were conserved locally (cf. Fig. 7, "decoupled").

Since $1 - \Gamma \not\ll 1$, the Keplerian angular velocity $\omega_{\rm Kep} = \sqrt{GM/R^3} \propto$

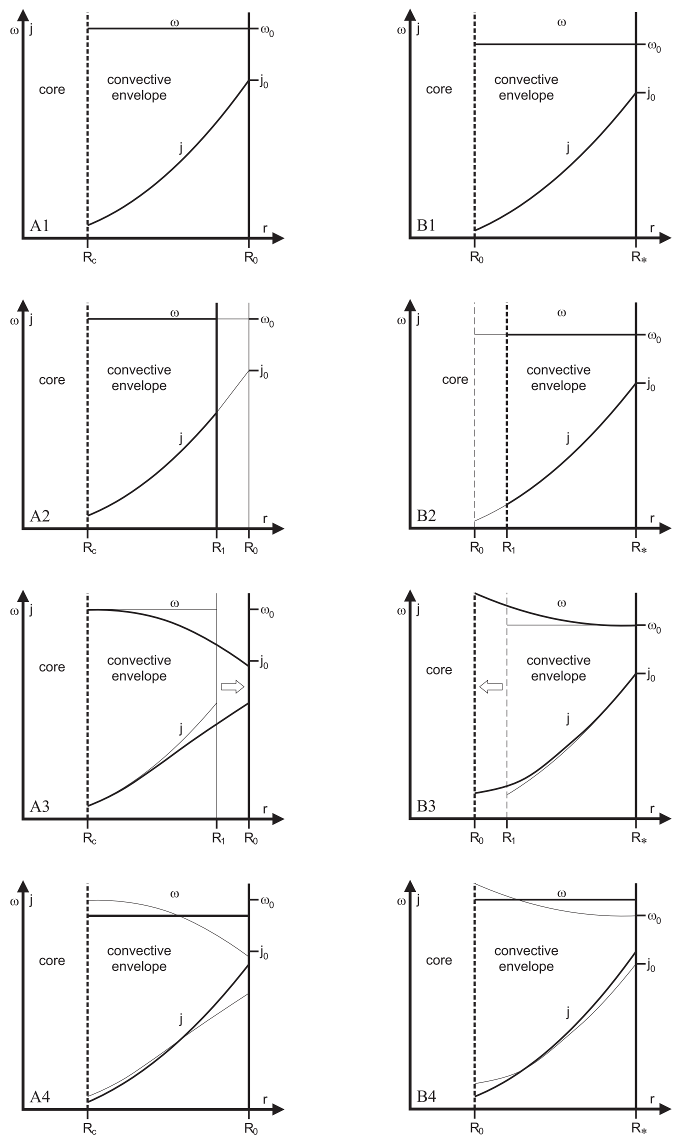

Figure 6. Mass loss from a rigidly rotating stellar envelope from the surface (case A: left panels) and through its lower boundary (case B: right panels). The continuous process is split up into three steps. First (panels 1 → 2), mass gets lost from the envelope, secondly, the envelope restores its original (inner or outer) radius R_0 (panels 2 → 3) by expansion, and third (panels 3 → 4) the specific angular momentum j is redistributed such that rigid rotation (i.e. $\omega(r)$ =const.) is restored. This leads to spin-down (spin-up) and decrease (increase) of the mean specific angular momentum for the case of mass loss through the upper (lower) boundary of the rigidly rotating stellar envelope. Thin lines show the state of the preceding step.

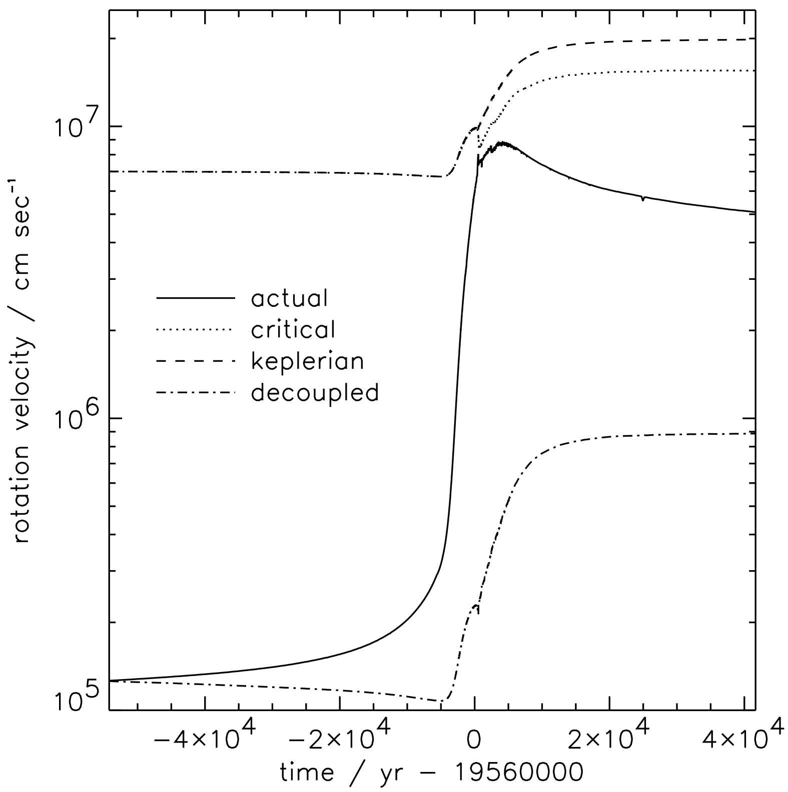

Figure 7. Equatorial rotation velocity as a function of time (solid line) compared to the Keplerian (dashed line) and the critical (dotted line) rotation rate; the latter two are different by the factor $1-\Gamma$. During the red supergiant phase it is $\Gamma \ll 1$ and the two lines coincide, while during the blue supergiant phase Γ rises to 0.4. The dash-dotted line shows the evolution of the surface rotation rate if there were no angular momentum transport in the convective envelope.

$f^{3/2}$ can be used as a rough approximation for the critical rotation frequency $\omega_{\rm c}$. Thus, Ω scales as $f^{1/2}$, and the mere contraction, without any angular momentum transport, would bring the star by a factor ~ 3 closer to its critical rotation rate. The numerical value of $\Gamma \approx 2\,10^{-3}$ on the RSG found in our calculation and $\Gamma \approx 0.4$ on the BSG would lead to an increase of Ω by ~ 3.9 from the RSG to the BSG, again assuming that j were conserved locally. Actually, the star gets from $\Omega \approx 0.01$ on the red supergiant branch at the beginning of central helium burning to critical rotation ($\Omega = 1$) during the RSG/BSG transition, even though a considerable loss of angular momentum due to stellar wind mass loss is included in our model (cf.

Fig. 7).

In fact, the model sequence would have exceeded the Ω-limit if we would not have applied a mass loss increase for $\Omega \rightarrow 1$ as in the massive main sequence models described in Section 2. I.e., like in that case, we find that stars undergoing a blue loop are likely to arrive at critical rotation, and they may consequently develop a slow equatorial outflow, which makes them candidates for B[e] stars. However, in contrast to the very massive star models discussed in Sect. 2, the B[e]-phase according to the blue loop scenario is much shorter, i.e. only some 10^4 yr, and correspondingly the amount of mass lost during that phase is much smaller. We want to mention here at least one star of which we know that it performed a red-blue evolution and for which the effect described here almost certainly played a role: the progenitor of Supernova 1987A (cf. Arnett et al. 1989, Langer et al. 1989; cf. also Woosley et al. 1997), and the possible twin of it as described by Brandner et al. (1997).

4. R4 and the binary scenario

In this Section, we want to investigate a binary scenario as possible explanation of the B[e]-phenomenon, which has emerged while trying to understand the properties of the B[e] supergiant R4, on which Franz-Josef Zickgraf has drawn our attention, and which is investigated in detail by Zickgraf et al. (1996b). These authors find that R4 is a spectroscopic binary with a period of 21.3 yr and an orbital separation of 23 A.U. ($\sim 5000\,R_\odot$). Presently, the system consists of a B[e] star with a luminosity of $L = 10^5\,L_\odot$ and an evolved ($T_{\rm eff} \simeq 9500\,$K) A star with $L = 1.4\,10^4\,L_\odot$.

These parameters lead immediately to an apparent contradiction: since the A star has already evolved off the main sequence, the almost 10 times more luminous companion should have become a supernova long time ago. Also, the orbital separation is too large in order to solve this contradiction by assuming that the A star was once more massive and shed mass onto the B component.

However, another binary scenario may be able to not only explain the properties of both stars but also the existence of a disk around the B star. In this scenario, the system was initially a triple system consisting of a close pair of stars with an initial mass of $\sim 10\,M_\odot$ for each of them, and another $\sim 10\,M_\odot$ star surrounding the pair on a wide orbit. The latter star is identical to the present A star, and only serves as a clock without ever interacting with the close pair.

The evolution of the close pair is sketched in Fig. 8. In order to end the evolution with the merging of both stars, the more massive component — the primary — must lose mass to the secondary at a high rate by Roche-

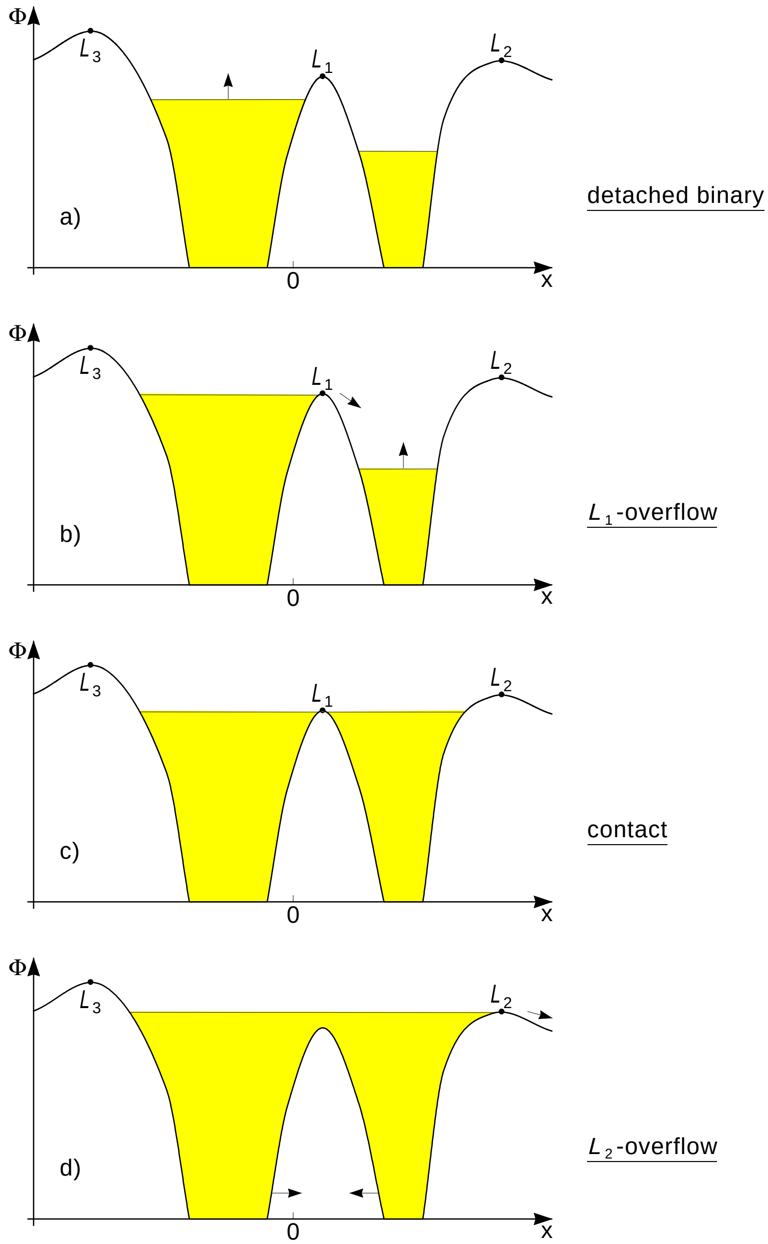

Figure 8. Schematic evolution of the spatial extents of two stars of a close binary system which evolves through a contact phase and develops mass outflow through the second Lagrange-point ("L_2-overflow"), which can be responsible for the formation of a circumsystem disk or ring (see text for details).

lobe overflow through the first Lagrange-point L_1. If the accretion time scale $\tau_{\dot{\rm M}} = M/\dot{M}$ of the secondary is shorter than its thermal time scale $\tau_{\rm KH} = GM^2/(RL)$ it will swell and start filling its own Roche-lobe. As the expansion will not stop here, matter will soon leave the system through the second Lagrange-point L_2, and the corresponding angular momentum loss will bring both stars continuously closer, with the merging of both stars as final result. Since the outflowing matter leaves the system in the orbital plane, i.e. through the L_2-point which rotates around the center of mass, it may form a ring or disk-like structure.

The most favorable initial conditions for such a scenario may be that the mass transfer process starts after the primary exhausted hydrogen in its center, but shortly before its subsequent expansion leads to the formation of a convective envelope (i.e., late Case B). Would the mass transfer start much earlier, i.e. during the core hydrogen burning phase of the primary (Case A), the merger star, which is now much more massive than the distant companion, would evolve into a supernova before the latter can leave the main sequence. Would the mass transfer occur when the primary has a red supergiant structure (Case C), both stars might merge as well, but so quickly that a common envelope is formed which, if at all, might tend to leave the star rather in a spherically symmetric way.

The merger remnant of a Case B binary would in fact, due to its small helium core mass, evolve into a hot blue supergiant star close to the main sequence in the HR diagram (cf. Podsiadlowski et al. 1992, Braun & Langer 1995), which fits to the effective temperature of the B[e] component of R4 of 27 000 K. Due to a strong helium overabundance in its envelope, it would appear overluminous for its mass, which seems actually to be observed: Zickgraf et al. (1996b) find a present mass of $\sim 13\,M_\odot$, but derive an initial mass of $\sim 20\,M_\odot$ from the comparison of its luminosity with standard single star evolutionary tracks.

5. Conclusions

In the previous Sections, we have presented three different evolutionary scenarios for the formation of a disk-like structure around a hot evolved massive star. The properties of the corresponding B[e] candidate models are compared in Table 1.

First, we can compare the positions of our B[e] candidates with the observed distribution of Magellanic Cloud B[e] supergiants in the HR diagram (Gummersbach et al. 1995; cf. also de Winter & van den Ancker 1997, and Zickgraf, this volume). Clearly, our main sequence scenario (Sect. 2) would not work for the stars with luminosities below $\sim 10^5\,L_\odot$. However, those stars are found to be comparatively cool ($\log T_{\rm eff} \simeq 4.1$) so that they fit

TABLE 1. Comparison of observable properties predicted by the three B[e] evolutionary scenarios discussed in this paper.

	very massive main sequence star at the Ω-limit	supergiant on blueward excursion from Hayashi line	single star remnant of binary merger
$T_{\rm eff}$	ok for most luminous B[e]s	ok for less luminous B[e]s	scatter in $T_{\rm eff}$ plausible
luminosity	ok for most luminous B[e]s	ok for less luminous B[e]s	scatter in L plausible
time scale	some 10^5yr	some 10^4yr	(?)
time integr. disk mass	$\sim 5\,M_\odot$	$\sim 0.1\,M_\odot$	$\sim 5\,M_\odot$(?)

quite well to the blue loop scenario outlined in Sect. 3. The very luminous group of B[e] stars investigated by Gummersbach et al., which can not correspond to stars on a blue loop, are mostly rather hot ($\log T_{\rm eff} \simeq 4.4$) and might well be main sequence stars. I.e., both scenarios together cover roughly the range of observed B[e] supergiants in the HR diagram.

For the binary scenario (Sect. 4), there are no quantitative models, but it appears likely that, assuming a scatter in the initial orbital parameters and in the initial mass ratio, the observed scatter in the HR diagram could also be reproduced.

As indicated in Table 1, the time scale of the B[e] phenomenon and the time integrated equatorial mass loss are rather different for the three models. While the main sequence model predicts a long B[e] life time with several solar masses expelled through the disk wind, the blue loop scenario predicts a shorter life time and less equatorial mass loss. The expected number of B[e] stars from both scenarios may still be comparable due to the steep decline of the initial mass function for larger masses.

B[e] time scale and disk mass loss are least clear for the binary scenario. During the L_2-overflow phase (cf. Fig. 8) certainly of the order of several solar masses of matter are lost. However, it is unclear which fraction of that is pushed to infinity and which fraction forms a disk. Another possibility which might be considered is that the star which is formed by the merger process is an extremely rapid rotator, and consequently has an equatorially focussed wind. In this case, the time scale of the B[e] phenomenon would be of the order of the spin-down time scale of the star.

Finally, we may also consider to combine the three scenarios discussed above. For example, if the merger of two stars is a red supergiant, e.g. in a Case C system, it may have considerably more angular momentum than red supergiant formed from single stars. If such a star evolves into a blue supergiant, our blue loop scenario must operate again, with an even stronger effect than in the single star case worked out in Section 3. It may be interesting to note in this context that Podsiadlowski (1997) suggested such a merger scenario for the presupernova evolution of SN 1987A.

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